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THERMAL EVOLUTION OF ISOLATED NEUTRON STARS

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Abstract. I give a general overview of the theory of neutron star cooling, emphasizing the intuitive understanding of the effects of the various physical ingredients, including the recently proposed accreted envelopes and neutrino emission from Cooper pair breaking and formation. I describe how the present data may be compatible with both the 'standard model' and the fast cooling scenarios, with or without invoking extensive presence of baryon pairing.

1. Introduction

Understanding the interior of neutron stars is a challenge to human intelligence and the study of their thermal evolution is one of the few possible methods through which such understanding (or misunderstanding) can be confronted with observations. The recent flow of data from *ROSAT* and the expected future avalanche of high quality data from *XMM* and *AXAF*, among others, has given a strong impetus to the field, which has resulted in many new important developments. My purpose here is to present the basic picture, emphasizing an intuitive understanding of the effect of the most important physical ingredients involved in this study. I have tried to refer to most of the recent works in the field in order to provide the reader with key entries into the literature.

2. Basic Physics of Neutron Star Cooling

The thermal evolution of a neutron star is basically given by the energy conservation equation

$$\frac{dE_{th}}{dt} = -L \qquad (1)$$

where E_{th} is the thermal energy content of the star and L the total luminosity, supplemented by the heat transport equation (and general relativistic corrections) [44]. L naturally divides into the surface photon emission $L_\gamma = 4\pi R^2 \sigma T_e^4$ and the internal neutrino emission L_ν. Internal heating mechanisms can be included in L as negative energy sinks H:

$$L = L_\gamma + L_\nu - H. \tag{2}$$

The following subsections give a brief description of the most important physics ingredients needed, whose recipes are put into an appropriate stellar evolution code [25, 24] which is easily crunched by a workstation.

2.1. THE EQUATION OF STATE

The internal structure of the neutron star is determined by the equation of state (EOS) which gives us two different inputs: the radial density profile and the 'chemical' composition, i.e., the type of particles present. I will use the classical Friedman & Pandharipande (FP) EOS [6] to determine the density profile and will add the core structure 'by hand': in principle, this is not the best thing to do, but there is so much uncertainty regarding the other ingredients described below that using EOSs appropriate to each scenario would not make much difference for this general description of neutron star cooling, and the difference could always be canceled by slightly adjusting another ingredient.

2.2. THE ENVELOPE

Traditionally considered to be the layer beneath the atmosphere down to a boundary density of $\rho_b = 10^{10}$ gm cm^{-3}, the envelope is a region of utmost importance. Its temperature gradient is very large and when the interior is isothermal it is the only region where a temperature gradient still persists. The envelope has generally been considered to be formed of iron-like nuclei and in this case, as a rule of thumb, the temperature at its base T_b and the effective temperature T_e are related by [10]

$$T_e \overset{\sim}{\propto} T_b^{0.5} \qquad\qquad T_e \approx 10^6 \text{ K} \longleftrightarrow T_b \approx 10^8 \text{ K}. \tag{3}$$

When the star is isothermal T_b is equal to the interior temperature.

Recently, Chabrier, Potekhin & Yakovlev [4] (see also [30, 39]) showed that the presence of light elements in the envelope strongly affects the heat transport and results, for a given T_b and at not too low a temperature, in higher surface temperatures compared to the previous models with beta-equilibrium matter. *This strongly affects the cooling and can change radically the conclusions drawn by comparing models to the data.*

The pulsar magnetic field, if higher than about 10^{11} G, also seriously affects the envelope, but, when realistic surface field configurations are considered [26, 38], the overall effect is not as large as has sometimes been claimed, at least in the case of iron envelopes; the case of magnetized accreted envelopes remains to be studied.

2.3. NEUTRINO EMISSION

The dominant neutrino emission processes occur in the core. If neutrons and protons are the only baryons present and the proton fraction is not too high, we are within the 'standard model' with the modified Urca process and its associated bremstrahlung brothers giving a neutrino emissivity

$$\epsilon_\nu^{MU} \;\cong\; 10^{20-21} \cdot T_9^8 \;\; \mathrm{erg\ cm^{-3}\ s^{-1}}. \tag{4}$$

Any change to this standard picture will almost always increase the neutrino emission by orders of magnitude: a larger proton fraction (which makes the direct Urca process possible [13]), a pion or kaon condensate, hyperons, quarks, etc... (see [29, 31] for reviews). It is thus convenient to divide the many possible scenarios into *slow* and *fast* neutrino cooling scenarios [21], the former being the 'standard model' and the latter being any of the other ones [1]. I will summarize the emissivity of the fast neutrino emission processes as

$$\epsilon_\nu^{\mathcal{N}} = 10^{\mathcal{N}} \cdot T_9^6 \;\; \mathrm{erg\ cm^{-3}\ s^{-1}} \tag{5}$$

with $\mathcal{N}$ ranging from about 24 (kaon condensate) up to 27 (direct Urca) [21] and a T^6 dependence instead of T^8 as in Eq. 4.

Besides the controversy about which - if any - fast neutrino emission process is allowed, there is still considerable uncertainty on the modified Urca rate. Voskresensky & Senatorov [50] proposed that medium effects may increase it by up to $2-3$ orders of magnitude, a possibility which obviously has a strong impact on the cooling [36].

Although proposed many years ago [5, 51], the neutrino emission by neutron (and proton) Cooper pair breaking and formation (PBF) has mysteriously been completely neglected until very recently [36]. Its neutrino emissivity can be comparable to or even higher than the modified Urca value but it only acts at $T < T_c$ and is rapidly suppressed when $T \ll T_c$: the result is a pulse of neutrino emission and enhanced cooling of the layer undergoing the pairing phase transition until $T \ll T_c$.

Neutrinos are also emitted in the crust, but their effect is practically negligible except at the early stages (~ 100 yrs).

[1] I like to distinguish *scenario* from *model*, the latter being a particular realization of the former which the computer can crank with all its details. Nevertheless, I will still use the term the 'standard model' instead of 'standard scenario'.

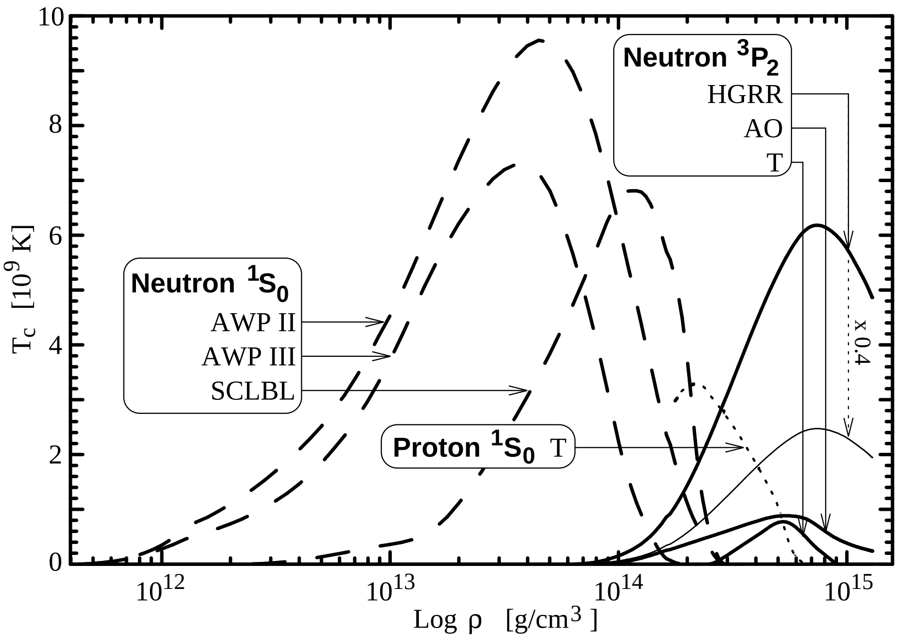

Figure 1. **Pairing critical temperatures used in this paper.** Neutron $^1S_0$: AWP II & AWP III from [1], and SCLBL from [37] which are the most reliable values calculated to date. Neutron $^3P_2$: HGRR from [11], and the same scaled down by 0.4, AO from [3], and T from [41]; the last two being more realistic than the first one. Proton $^1S_0$: T from [42] which is close to the better calculation of [52]. The maximum density shown is the central density of my 1.4 $M_\odot$ FP neutron star.

2.4. SUPERFLUIDITY AND SUPERCONDUCTIVITY

In the pairing of the neutrons (= superfluidity) and protons (= superconductivity), the Cooper pairs form in some angular momentum state which is usually $^1S_0$ at low momentum and a mixed $^3P_2 - {}^3F_2$ state at higher momentum [33]. As shown in Figure 1, there are still large uncertainties on the values of the critical temperatures T_c at which these pairing phase transitions occur, particularly in the core. The pairing phenomenon introduces a gap Δ in the single particle excitation spectrum which, when appropriately defined, is related to T_c by $\Delta(0) \simeq 1.75 T_c$. At $T \ll T_c$ this gap produces a strong suppression of the specific heat of the paired component and of the neutrino emission processes in which this component participates:

$$c_v^{(N)} \;\to\; c_v^{(P)} = \mathcal{R}_c \cdot c_v^{(N)} \qquad\qquad \epsilon_\nu^{(N)} \;\to\; \epsilon_\nu^{(P)} = \mathcal{R}_\nu \cdot \epsilon_\nu^{(N)} \tag{6}$$

where 'N' stands for 'normal' and 'P' for 'paired'. The suppression coefficients $\mathcal{R}_c$ and $\mathcal{R}_\nu$, have been described in detail by Levenfish & Yakovlev [15, 16, 53].

2.5. INTERNAL HEATING

Several mechanisms of internal heating have been proposed, for example: friction due to the differential rotation of the crustal neutron $^1S_0$ superfluid [2]; dissipative processes due to the core proton $^1S_0$ superconductor vortex lines [34]; release of 'chemical' energy due to the readjustment of the chemical equilibrium of the core induced by the spin-down of the pulsar [32]. The first of these is potentially the most efficient for young pulsars and has already been studied in some detail [40, 46, 48]. I will adopt it in a simple version [23], writing the heating rate as

$$H(t) = J_{44} \cdot 10^{40} \cdot \left(\frac{t + \tau_0}{100 \text{ yrs}} \right)^{-3/2} \quad \text{erg s}^{-1} \tag{7}$$

where t is the pulsar age, $\tau_0 = 300$ yrs a typical spin-down time scale, and J_{44} the differential angular momentum of the frictionally coupled crustal neutron superfluid in units of 10^{44} g cm^2 rad s^{-1}. This expression assumes a standard spin-down rate from magnetic dipole radiation and is similar to the one used in [40, 46]. This heating is distributed within the superfluid layers of the inner crust. I adopt the value $J_{44} = 3.1$ which corresponds to moderately strong heating, compatible with the size of the crust for the FP EOS. Very different heating rates are of course possible [48].

3. An overview of the various scenarios

As the previous section made clear, we have a plethora of effects which may affect the thermal evolution of a neutron star. Several of them have been studied in detail in the literature while others have been proposed, or simply unearthed, very recently and need more consideration. I will show here a series of cooling curves which illustrate these various effects in order to see them at work and try to understand them intuitively. Figures 2 to 4 should be deciphered by careful comparison with Figure 1. Fast neutrino emission, when used, is plugged in at densities $\rho > \rho_{cr} = 1.1 \cdot 10^{15}$ gm cm^{-3}.

3.1. THE GENERAL PRE-1996 PICTURE

I show in Figure 2 a set of curves to illustrate the main cooling scenarios and the effect of baryon pairing. At an early stage ($t \lesssim 10 - 100$ yrs) all models have the same surface temperature: the surface does not know yet what is happening in the core, because of the finite time scale for heat diffusion through the crust [19], and its temperature is controlled by plasma neutrino emission in the outer crust and by surface photon emission. All models use the same FP EOS and thus have the same crust and hence exactly the same surface temperature at this stage.

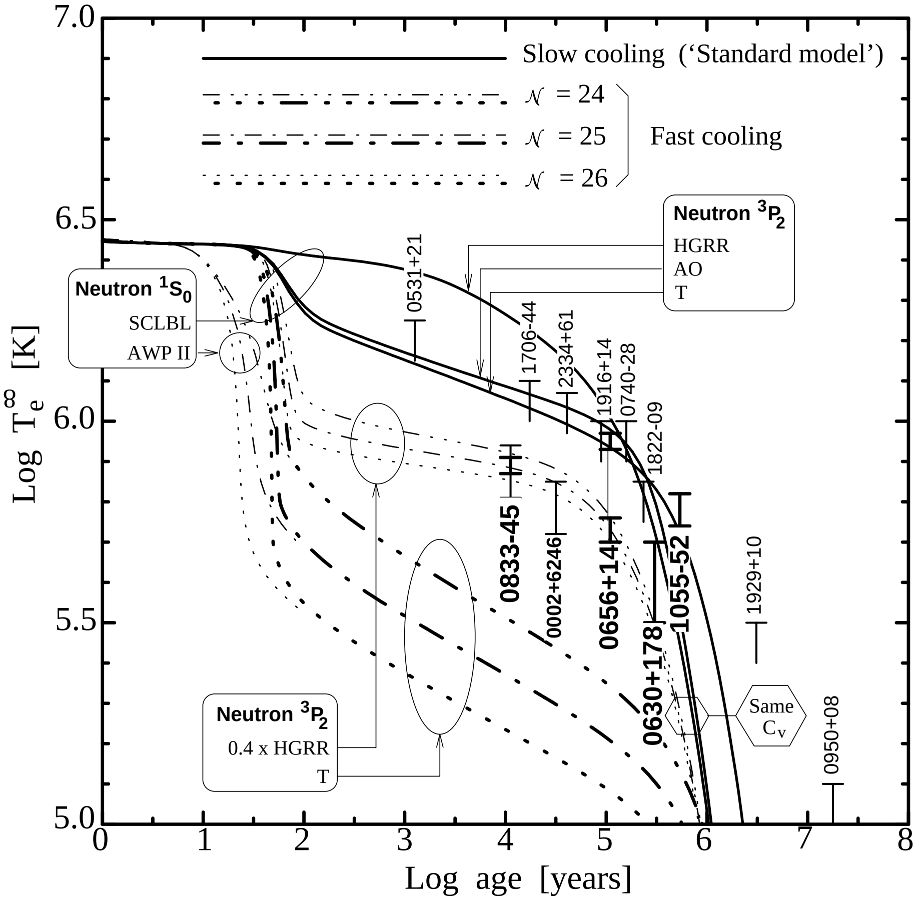

Figure 2. **Typical behavior of slow ('standard') and fast cooling scenarios.** 1.4 $M_\odot$ neutron stars with the FP EOS. The cases with $\mathcal{N}$ = 24, 25, and 26 (see Eq. 5) correspond approximately to the effect of a kaon condensate, pion condensate, and the direct Urca process (with hyperons or nucleons), respectively. The various curves, within each scenario, show the effect of various assumptions about pairing, following the notations of Figure 1: all models use the proton $^1S_0$ T_c 'T', and the neutron $^1S_0$ and $^3P_2$ T_c's are as labeled. All models have non-magnetized iron envelopes from [4]. Neither PBF neutrino emission nor heating are included.
The main effect of pairing in the crust (neutron $^1S_0$) is to shorten the length of the early plateau. Core pairing suppresses the neutrino emission, which results in a higher T_e during the neutrino cooling era (age from $\sim$ 100 to $\sim 10^5$ yrs), and the specific heat, which results in faster cooling during the photon cooling era (age above $\sim 10^5$ yrs). The reduction of the specific heat during the neutrino cooling era does not show up as much as during the photon cooling era due to the small slope of the curves at this phase.
The references to the data can be found in [23]: in short, the bigger the label the better the data. All points are really upper limits (in several cases based on a non-detection of the pulsar) but for the radio pulsars 0833-45 (Vela), 0656+14, 0630+178 (Geminga) [12], 1055-52, and the neutron star 0002+6246, there is good evidence that the observed X-rays are from surface thermal emission. Uncertainty on the temperature estimate is illustrated in the case of PSR 0656+14 where two values are reported. For more details see [43, 23].

After this early *plateau* the surface temperature decreases more or less strongly, depending on the evolution of the core: this is the *isothermalization phase* during which the heat content of the crust (and outer core) flows into the core, where it is evaporated in neutrinos. At the end of this phase the temperature profile inside the star is flat, i.e., the star is isothermal, except for the strong temperature gradient within the envelope. The time needed for isothermalization is $\approx D^2 \cdot C_v/\lambda$ where D is the thickness of the insulating layer (= crust + outer core), C_v its specific heat, and λ its thermal conductivity. Observation of a neutron star at this stage would directly 'measure' the thickness D [14]. However, this requires accurate pairing gaps in the crust to calculate C_v since pairing reduces it. The two sets of fast cooling curves in Figure 2 with the neutron $^1\mathrm{S}_0$ pairing from AWP II and SCLBL show it clearly: in the AWP II case the gap extends to lower densities, implying a lower specific heat and thus an earlier temperature drop. Moreover, neutrino emission by the PBF process (see § 3.2.2) complicates the situation. In the case of a strange star [9], this plateau lasts only about a year [20] due to the absence of an inner crust, i.e., here D is very small.

Once the star is isothermal we can distinguish a *neutrino cooling era* followed by a *photon cooling era* depending on which energy loss mechanism is driving the evolution: the change from the former to the latter shows itself as a change in the slope of the cooling curves, occuring around $t \sim 10^5$ yrs.

During the neutrino cooling era we see a clear difference between the 'standard model' and the fast cooling scenarios. Within the fast cooling scenarios, the cases with the neutron $^3\mathrm{P}_2$ T_c from 'T' [41] show the fast cooling unsuppressed: this gap, as well as the proton gap, vanishes within the inner core where fast neutrino emission occurs (Figure 1). The resulting surface temperatures are much below the observed ones and it had been proposed [24, 25] that pairing gaps extending down to the center of the star could control the fast neutrino emission and keep the star much warmer. The suppression of the neutrino emission by the gap is of course very sensitive to T_c, higher T_c's resulting in higher surface temperatures [25, 24]. One could actually use this to 'measure' T_c in the core of a neutron star by fitting the cooling curves to the data [8, 17, 22, 24]. This is what I have done in the models with neutron $^3\mathrm{P}_2$ T_c labeled as '0.4 x HGRR': comparison with Figure 1 shows that T_c's around 2×10^9 K are needed. This 'measurement' needs exact suppression factors (Eq. 6) as used recently in [7, 8, 17, 22, 23] and implies values of T_c higher than when simple Boltzmann factors are used as, recently, in [24, 25, 35, 36, 45, 47]. Notice that when the fast neutrino emission is suppressed by pairing, the resulting cooling is only very weakly dependent on the exact cooling agent (here the value of $\mathcal{N}$), the EOS, and the critical density at which this agent starts operating [22]. Within the 'standard model' the effect of pairing is similar

but less spectacular: the curve with the neutron ${}^3\mathrm{P}_2$ T_c from 'AO' [3] is slightly warmer than the one with 'T' [41] since the latter T_c is lower and also vanishes in the inner core, while the curve with the T_c 'HGRR' [11] is much warmer because the neutrino emission has been suppressed very early on, in the whole core, due to the high value of T_c.

During the photon cooling era the temperature decreases much faster (on a log scale) and the different cooling scenarios largely overlap: observations of neutron stars of age $\gtrsim 10^5$ yrs do not allow us to distinguish between them [21]. The cooling rate now depends crucially on the total specific heat (and the structure of the envelope, as shown in § 3.2.1). For example, the two 'standard' cooling curves 'HGRR' and 'AO' merge at this stage, despite the very different T_c's, because in both cases the neutrons are paired in the whole core and $T << T_c$, while the 'T' curve cools more slowly since its inner core neutrons are not paired and its specific heat is thus larger. Moreover, notice that the fast cooling models with the '0.4 x HGRR' T_c asymptotically approach the 'standard' cooling curves 'HGRR' and 'AO': at this time they all have the neutrons paired in the whole core and thus the same specific heat.

It is also most instructive to look at detailed temperature profiles for the various scenarios: many have been published, for example in [14, 18, 19, 46, 49] and the models of this paper are available on the Web.

Magnetic fields in the envelope do alter the picture somewhat, but not strongly [26, 38]: light elements in the envelope and PBF neutrino emission have much more dramatic effects as shown in the next section.

3.2. NEW DEVELOPMENTS

The picture presented in the previous section may change dramatically due to three new ingredients introduced recently: the possible presence of light elements in the upper layers of the neutron star, the neutrino emission by the breaking and formation of Cooper pairs, and the possibility of substantial in-medium enhancement of the modified Urca process.

3.2.1. *Accreted envelopes (and heating)*

The light elements in the envelope increase the heat transport, resulting in a higher T_e for a given T_b, as long as T_e is not too low [4, 30, 39]. During the early stage and the neutrino cooling era the surface temperature simply follows the evolution of the interior temperature and does not affect the cooling rate. As a result, at these stages a neutron star with an accreted envelope will undergo an evolution parallel to the evolution of the same neutron star with an iron envelope but with a higher T_e. When photon cooling takes over the situation is reversed, since for a given T_b (i.e., in-

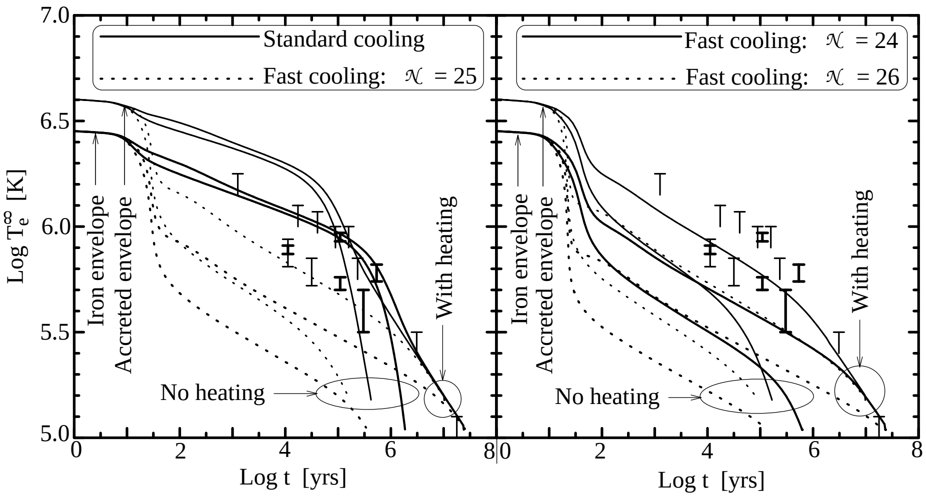

Figure 3. **Effect of an accreted envelope and heating on the cooling.** 1.4 $M_\odot$ neutron stars star with the FP EOS. The fast neutrino emission is plugged in as described in § 2.3. All models use T_c's from 'AWP II' [1] and 'T' [41] for neutrons, and 'T' [42] for protons (see Figure 1): in these models there is *no* pairing in the inner core. The presence of light elements in the envelope raises the surface temperature during the neutrino cooling era and then hastens the cooling during the photon cooling era. The internal heating works at all times but is of course more efficient when the heat content, i.e., interior temperature, of the star is low. When the initial heat content of the star has been completely lost the cooling is independent of the previous history and envelope structure since the luminosity is simply equal to the heating rate. Notice that the temperature estimate of Geminga had been argued to require extensive pairing in the core to be compatible with cooling models [21], both 'standard' and fast: in models with accreted envelopes, as shown here and in [4, 30], *extensive baryon pairing in the core is not necessary*. The data points are labeled in Figure 2.

terior temperature and thus specific heat) the surface emission is stronger in the case of an accreted envelope because T_e is higher. Thus, the cooling trajectories of these two neutron stars cross and the temperature of the one with an accreted envelope drops much faster. This behavior is illustrated in Figure 3 above and, even more clearly, in the Figure 2 of [39].

If, moreover, internal heating is included, one can obtain higher temperatures at all ages (nothing new in this of course, see [40, 46, 48]). This is extremely important for the fast cooling scenarios for which the results of Figure 3 show that *it is possible to obtain temperatures in agreement with the presently available data with models where the gap vanishes within the inner core* [23], in contradistinction to the results of § 3.1. The case $\mathcal{N}$ = 24 (i.e. a kaon condensate or a pion condensate with its neutrino emission reduced by medium effects [45]) appears to be the most favorable but even the direct Urca process ($\mathcal{N}$ = 26 - 27) could be acceptable.

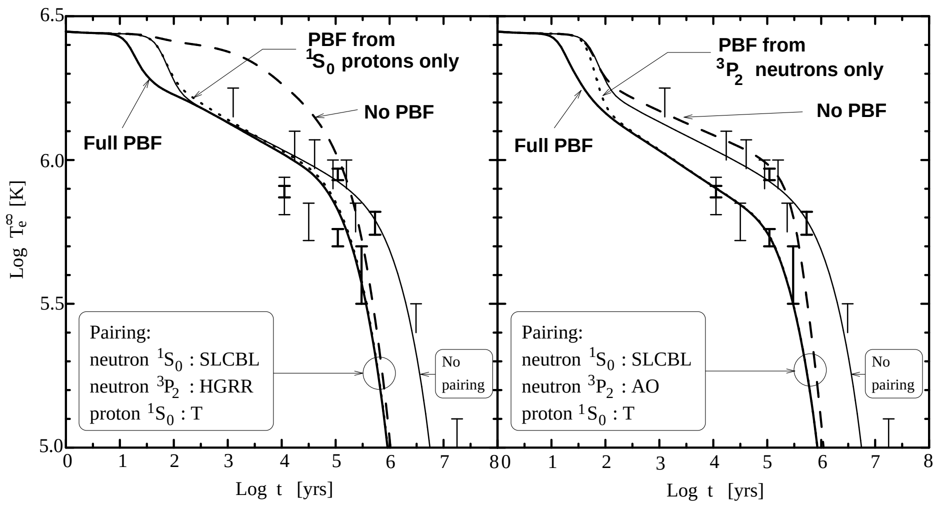

Figure 4. **Effect of the neutrino emission by the Pair Breaking and Formation (PBF) process on the 'standard' cooling.**
1.4 $M_\odot$ neutron stars with the FP EOS. The pairing T_c's are as labeled (see Figure 1). These models have an iron envelope and no heating is included.
The PBF emission in the crust hastens the isothermalization but has little effect later on. Once the star is isothermal, the cooling may be driven by the PBF emission from either core proton $^1S_0$ pairing (left panel) or core neutron $^3P_2$ pairing (right panel): it is the component with the lowest T_c which dominates since its PBF neutrino emission happens later. In these particular choices of pairing, the neutron $^3P_2$ PBF emission in the right panel is more efficient than the proton $^1S_0$ PBF emission in the left panel since a larger mass of paired matter is involved. The data points are labeled in Figure 2.

3.2.2. *Neutrino emission from Cooper pair breaking and formation*

Besides its suppressive effect, the occurrence of pairing does induce a strong, but short lived, neutrino emission [5, 51] due to the constant breaking and formation of Cooper pairs (§ 2.3). I show in Figure 4 'standard' cooling curves with these PBF mechanisms included: the results are quite sensitive to the values of T_c for neutrons and/or protons in the core but the effect can be sufficient to lower the star temperature significantly. One sees that, *within the 'standard' cooling scenario with the (also 'standard') PBF neutrino emission taken into account there may be no need of any 'exotic' process* (see Schaab *et al.* [36]). Notice, however, that none of the models shown in Figure 4 include any heating or an accreted envelope: inclusion of these could spoil the apparent agreement with the data.

3.2.3. *Enhanced modified Urca process*

If, as Voskresensky & Senatorov [50] proposed, the modified Urca rate is much more efficient than usually assumed, then all 'standard' cooling curves

are pushed down and *all data, as presently interpreted, could be compatible with the 'standard model'* (see Schaab *et al.* [36]). The actual efficiency of the modified Urca process is a very delicate, and controversial, problem but if neutron (or proton) pairing is present within the whole core it would be suppressed anyway and possibly still result in too high temperatures.

4. Conclusions

Apparently, our misunderstanding of the neutron star interior has progressed as much as our understanding. The present data seem to be compatible with both the standard scenario and the fast cooling scenarios, depending on the other assumptions made in building the models. However, the situation is far from desperate:

– On the observational side, future - and multiwavelength - observations will certainly be able to determine the chemical composition of the atmosphere [28], and thus probably of the envelope, of cooling neutron stars. Observations of old pulsars will refute or confirm the presence of internal heating [27] and help us to pin down its nature [48].

– On the theoretical side, what is badly needed, but yet can reasonably be expected to be obtained, is a modified Urca rate that everybody will agree upon (at least the order of magnitude) and reliable calculations of the pairing gaps for neutrons and protons in 'standard' matter. This will allow us to make more definite predictions for the standard scenario. Moreover, much work is constantly being done on all the fast cooling scenarios.

Finally, the good news is that, thanks to the accreted envelopes, the fast cooling scenarios are viable without having to invoke the doubtful presence of pairing at the extreme densities reached in 'exotic' neutron stars [23].

Data files for all cooling curves shown here are available on the Web at:
http://www.astroscu.unam.mx/neutrones/NS-Cooler/NS-Cooler.html

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References

1. Ainsworth, T.L., Wambach, J. & Pines, D. 1989, Phys. Lett., B222, 173
2. Alpar, M.A., Anderson, P.W., Pines, D., & Shaham, J. 1984, ApJ, 276, 325
3. Amundsen, L. & Østgaard, E. 1985, Nucl. Phys., A442, 163
4. Chabrier, G., Potekhin, A.Y., & Yakovlev, D. G. 1997, ApJ, 477, L99
5. Flowers, E., Ruderman, M., & Sutherland, P. 1976, ApJ, 205, 541
6. Friedman, B., & Pandharipande, V.R. 1981, Nucl. Phys., A361, 502
7. Gnedin, O.Yu., & Yakovlev, D.G. 1993, Astron. Lett., 19, 104
8. Gnedin, O.Yu., Yakovlev, D.G., & Shibanov, Yu.A. 1994, Astron. Lett., 20, 409
9. Glendenning, N. 1997, contribution to these proceedings

10. Gudmundsson, E.H., Pethick, C.J. & Epstein, R.I. 1983, ApJ, 272, 286
11. Hoffberg, M., Glassgold, A.E., Richardson, R.W. & Ruderman, M. 1970, Phys. Rev. Lett., 24, 775
12. Kuzmin, A.D., & Losovsky, B.Ya. 1997, Astron. Lett., in press (May 1997 issue)
13. Lattimer, J.M., Pethick, C. ., Prakash, M. & Haensel, P. 1991, Phys. Rev. Lett., 66, 2701
14. Lattimer, J.M., Van Riper, K.A., Prakash, M., & Prakash, M. 1994, ApJ, 425, 802
15. Levenfish, K.P., & Yakovlev, D.G. 1994a, Astron. Lett., 20, 54
16. Levenfish, K.P., & Yakovlev, D.G. 1994b, Astron. Rep., 71, 282
17. Levenfish, K.P., & Yakovlev, D.G. 1996, Astron. Lett., 22, 56
18. Muslimov, A., & Page, D. 1996, ApJ, 458, 347
19. Nomoto, K., & Tsuruta, S. 1987, ApJ, 312, 711
20. Page, D. 1992, in High Energy Phenomenology, eds M.A. Pérez & R. Huerta (Singapore: World Scientific) 347 (e-print: astro-ph/9602043)
21. Page, D. 1994, ApJ, 428, 250
22. Page, D. 1995, Rev. Mex. Fís., 41, Supl. 1, 178 (e-print: astro-ph/9501071)
23. Page, D. 1997, ApJ, 479, L43
24. Page, D. & Applegate, J. H. 1992, ApJ, 394, L17
25. Page, D. & Baron, E. 1990, ApJ, 354, L17; Erratum in ApJ, 382, L111
26. Page, D., & Sarmiento, A. 1996, ApJ, 473, 1067
27. Pavlov, G.G., Stringfellow, G.S., & Córdova, F.A. 1996, ApJ, 467, 370
28. Pavlov, G.G, Zavlin, V.E., Trümper, J., & Neuhäuser, R. 1996, ApJ, 472, L33
29. Pethick, C. J. 1992, Rev. Mod. Phys., 64, 1133
30. Potekhin, A. Y., Chabrier, G., & Yakovlev, D. G. 1997, A&A, in press
31. Prakash, M. 1994, Phys. Rep., 242, 297
32. Reisenegger, A. 1995, ApJ, 442, 749
33. Sauls, J. 1997, contribution to these proceedings
34. Sedrakian, A.D., & Sedrakian, D.M. 1993, ApJ, 413, 658
35. Schaab, Ch., Weber, F., Weigel, M.K., & Glendenning, N.K. 1996, Nucl. Phys, A605, 531
36. Schaab, Ch., Voskresensky, D., Sedrakian, A.D., Weber, F. & Weigel, M. K. 1997, A&A, in press
37. Schulze, H.-J., Cugnon, J., Lejeune, A., Baldo, M., & Lombardo, U. 1996, Phys. Lett., B375, 1
38. Shibanov, Yu. A., & Yakovlev, D. G. 1996, A&A, 309, 171
39. Shibanov, Yu. A., Potekhin, A.Y., Yakovlev, D.G., & Zavlin, V.E. 1997, contribution to these proceedings
40. Shibazaki, N., & Lamb, F.K. 1989, ApJ, 346, 808
41. Takatsuka, T. 1972, Prog. Theor. Phys., 48, 1517
42. Takatsuka, T. 1973, Prog. Theor. Phys., 50, 1754
43. Trümper, J. 1997, contribution to these proceedings
44. Tsuruta, S. 1986, Comments Astrophys., 11, 151
45. Umeda, H., Nomoto, K., Tsuruta, S., Muto, T., & Tatsumi, T. 1994, ApJ, 431, 309
46. Umeda, H., Shibazaki, N., Nomoto, K., & Tsuruta, S. 1993, ApJ, 408, 186
47. Umeda, H., Tsuruta, S., & Nomoto, K. 1994, ApJ, 433, 256
48. Van Riper, K. A., Link, B., & Epstein, R. I. 1995, ApJ, 448, 294
49. Van Riper, K. A. 1991, ApJS, 75, 449
50. Voskresensky, D. N., & Senatorov, A. V. 1986, Sov. Phys. JETP, 63, 885
51. Voskresensky, D. N., & Senatorov, A. V. 1987, Sov. J. Nucl. Phys., 45, 411
52. Wambach, J., Ainsworth, T.L., & Pines, D. 1991, in Neutron Star: Theory and Observation, eds. J. Ventura & D. Pines (Dordrecht: Kluwer Academic Publishers) 37
53. Yakovlev, D.G., & Levenfish, K.P. 1995, A&A, 297, 717

NOTE: four references not yet published are available on the Web:

[**9**] Glendenning, N. 1997, contribution to these proceedings
e-print: astro-ph/9706236

[**30**] Potekhin, A. Y., Chabrier, G., & Yakovlev, D. G. 1997, A&A, in press
e-print: astro-ph/9706148

[**36**] Schaab, Ch., Voskresensky, D., Sedrakian, A.D., Weber, F. & Weigel, M. K. 1997, A&A, in press
e-print: astro-ph/9605188

[**39**] Shibanov, Yu. A., Potekhin, A.Y., Yakovlev, D.G., & Zavlin, V.E. 1997, contribution to these proceedings.
Preprint at: http://stella.ioffe.rssi.ru/NSG/NSG-Pub1.html